THEORETICAL LINE PROFILES IN THE ULTRAVIOLET SPECTRA OF EARLY-TYPE STARS

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ABSTRACT

Theoretical ultraviolet spectra of two early B stars have been computed in detail from 911.6 to 3000 Å, including all lines expected to be wider than about 2 Å. The lines were assumed to be formed in pure absorption. Perturbations of both ions and electrons were included in the broadening of the hydrogen lines, while a Doppler-damping profile with a damping constant ten times the classical value was used for all other elements. Most of the lines are found shortward of 1410 Å where they have deep cores and many have extensive wings. These overlapping lines cause significant blanketing, amounting to at least 31 per cent of the total flux at $T_e = 25673°$ K and at least 20 per cent at $T_e = 31023°$ K. This effect must be included in the constant-flux criterion when computing the model atmospheres.

I. INTRODUCTION

Rocket-borne instruments soon will provide ultraviolet stellar spectra with resolutions of about 1 Å in which the strongest lines should appear. These spectra may contain many unexpected features, but nevertheless it is useful to make preliminary calculations of the lines with the best theory now available. The first rockets probably will be directed toward O and B stars, since the anticipated strong ultraviolet radiation from such hot atmospheres makes these types the most interesting and the easiest to observe.

Ultraviolet spectra of the Sun (Tousey 1963) have shown emission lines shortward of 1850 Å arising from the chromosphere and corona. It is not unreasonable that a hotter star also has a temperature inversion and outer layers where local thermodynamic equilibrium does not exist. However, in the Sun no chromospheric features are seen in the visible spectrum of the disk except in the cores of the very strongest lines, so that in an O or B star it would be surprising if non-equilibrium processes could enhance many lines above the ultraviolet continuum, which carries the bulk of the photon flux. For this reason all lines in this paper are assumed to be in absorption.

Gaustad and Spitzer (1961) have calculated the equivalent widths of the stronger ultraviolet lines to be expected in a uniform atmosphere with a temperature of 2×10^4 ° K and an electron pressure of 2×10^3 dyne cm^{-2}. Using the classical value for the damping constant they found forty-two lines wider than 1 Å and ten wider than 10 Å. With this constant ten times larger they estimated that the lines absorb 68 per cent of the flux below 1400 Å. This extreme blanketing effect indicates the importance of extending these calculations to obtain the blended line profiles in an atmosphere with a realistic temperature-pressure distribution. This paper reports the results of such a program.

Two models were chosen from a series of B-star model atmospheres by Miss Underhill (1962*a*). She used a surface gravity of 10^4 cm sec^{-2} and abundances by mass of 0.68 hydrogen and 0.32 helium, typical for main-sequence B stars. Electron scattering and the bound-free absorption of H I, H^-, He I, and He II were included in the opacity, but no line absorption. Table 1 lists the spectral types, boundary temperatures T_0, effective temperatures T_e, and electron pressures P_e at which $T = T_e$ for the two models. The spectral types were estimated from the calculated Balmer discontinuities. Strömgren (1964) compared the cooler model with his observed color index $u - b$, corresponding to a ratio of fluxes at 3500 and 4670 Å, and concluded that the spectral type is closer to B1 V. The hotter model also may be somewhat earlier in type. Recent models by Mihalas (1964) show excellent agreement with the temperature-pressure distributions adopted here except that the boundary temperatures are 1000° K lower. In the outermost layers, where the temperature drops rapidly due to the opacity in the Lyman continuum, Miss Underhill has very few integration points down to unit optical depth at wavelengths shortward of 911.6 Å.

II. THE LINE DATA

Table 2 contains the abundances by weight of the elements which are expected to have strong ultraviolet lines in these atmospheres. Helium and oxygen are added for reference, although probably none of their lines are wider than 2 Å. The abundances relative to oxygen are from the solar investigation of Goldberg, Müller, and Aller (1960) except for chlorine and argon which were obtained from studies of hot stars and listed by Aller (1958). These values were related to hydrogen by the solar oxygen-hydrogen ratio derived by Osterbrock and Rogerson (1961). A rediscussion of this ratio by Gaustad (1964) suggests that it and all the abundances except H and He in Table 2 might be reduced by a factor of $\frac{2}{3}$. The chlorine abundance determined from planetary nebulae is an additional factor of 5 lower than the one quoted here.

TABLE 1

CHARACTERISTICS OF THE MODELS

Spectrum	T_0 (° K)	T_e (° K)	P_e at $T=T_e$ (dyne cm^{-2})
B2 V	16000	25673	2.32×10^3
B0 V	21294	31023	2.14×10^3

TABLE 2

ABUNDANCES OF THE ELEMENTS BY WEIGHT

Element	Abundance	Element	Abundance
H	0.68	Si	8.8×10^{-4}
He	.29	S	6.7×10^{-4}
C	.0067	Cl	8.8×10^{-4}
N	.0014	Ar	6.3×10^{-4}
O	.015	Fe	2.2×10^{-4}
Al	0.45×10^{-4}		

The populations of the ion states of the elements were computed by Upson (1963) for the assumed atmospheres. Then the *Ultraviolet Multiplet Tables* of Moore (1950, 1952) were searched for the strongest lines of the abundant ions. Reference to the widths of Gaustad and Spitzer (1961) or to actual profile calculations showed that many of these lines were narrower than 1 Å and consequently were eliminated from this program which is aimed at studying only the most important ones. Table 3 contains those finally included in the blended profiles; it is hoped that for lines with available f-values none with total width at half-depth greater than 2 Å in these models has been omitted.

However, if such lines as N IV at 923.2 and Si IV at 1128.3 Å have f-values of the order of unity, they would be this wide. Other abundant ions which have strong ultraviolet lines in a list by Kelly are O IV, F III, Ne II, Ne III, Na III, Mg III, Al IV, Ar III, Fe IV, and Fe V. These should be considered when the multiplet designations and f-values are known.

In Table 3 the wavelengths and term designations are from the *Ultraviolet Multiplet Tables* of Moore (1950, 1952) while the excitation potentials are from her tables of

TABLE 3

THE STRONGEST ULTRAVIOLET LINES IN EARLY B STARS

Ion	Wavelength (Å)	Multiplet	J_l–J_u	Excitation (cm^{-1})	gf
L18	914.576	$1s\ ^2S$–$18p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.000540
L17	914.919	$1s\ ^2S$–$17p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.000642
L16	915.329	$1s\ ^2S$–$16p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.000770
L15	915.824	$1s\ ^2S$–$15p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.000938
L14	916.429	$1s\ ^2S$–$14p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.001154
L13	917.181	$1s\ ^2S$–$13p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.001446
L12	918.129	$1s\ ^2S$–$12p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.001842
L11	919.351	$1s\ ^2S$–$11\mathrm{p}\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.002402
L10	920.963	$1s\ ^2S$–$10p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.003210
L9	923.150	$1s\ ^2S$–$9p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.004432
L8	926.226	$1s\ ^2S$–$8p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.006366
L7	930.748	$1s\ ^2S$–$7p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.009628
L6 ϵ	937.804	$1s\ ^2S$–$6p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.01560
L5 δ	949.743	$1s\ ^2S$–$5p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.02788
L4 γ	972.537	$1s\ ^2S$–$4p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.05798
L3 β	1025.722	$1s\ ^2S$–$3p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.1582
L2 α	1215.670	$1s\ ^2S$–$2p\ ^2P^0$	$\frac{1}{2}$–$\frac{1}{2}$, $\frac{3}{2}$	0	0.8324
N II	915.603	$2p^2\ ^3P$–$2p^3\ ^3P^0$	0–1	0	0.47
	916.0	....	1–0, 1, 2	49.1	1.41
	916.700	....	2–1, 2	131.3	2.35
Ar II	919.78	$3p^5\ ^2P^0$–$3p^6\ ^2S$	$\frac{3}{2}$–$\frac{1}{2}$	0	1.72
	932.046	....	$\frac{1}{2}$–$\frac{1}{2}$	1432.0	0.86
C III	977.026	$2s^2\ ^1S$–$2p\ ^1P^0$	0–1	0	1.20
Cl IV	973.212	$3p^2\ ^3P$–$3p^3\ ^3D^0$	0–1	0	0.55
	977.7	....	1–1, 2	491	1.65
	984.952	....	2–3	1341	2.32
	985.749	....	2–1, 2	1341	0.44
N III	989.790	$2p\ ^2P^0$–$2p^2\ ^2D$	$\frac{1}{2}$–$\frac{3}{2}$	0	0.94
	991.514	....	$\frac{3}{2}$–$\frac{3}{2}$	174.5	0.19
	991.579	....	$\frac{3}{2}$–$\frac{5}{2}$	174.5	1.69
Si III	993.54	$3p\ ^3P^0$–$4s\ ^3S$	0–1	52630	0.20
	994.82	....	1–1	52758	0.60
	997.40	....	2–1	53019	1.00
Cl III	1005.280	$3p^3\ ^4S^0$–$3p^4\ ^4P$	$\frac{3}{2}$–$\frac{1}{2}$	0	0.56
	1008.777	....	$\frac{3}{2}$–$\frac{3}{2}$	0	1.10
	1015.023	....	$\frac{3}{2}$–$\frac{5}{2}$	0	1.66
C II	1009.854	$2p^2\ ^4P$–$2p^3\ ^4S^0$	$\frac{1}{2}$–$\frac{3}{2}$	43000.2	1.00
	1010.074	....	$\frac{3}{2}$–$\frac{3}{2}$	43021.8	2.00
	1010.369	....	$\frac{5}{2}$–$\frac{3}{2}$	43050.7	3.00
S III	1012.49	$3p^2\ ^3P$–$3p^3\ ^3P^0$	0–1	0	0.30
	1015.51	....	1–0, 1	297.2	0.52
	1015.76	....	1–2	297.2	0.38
	1021.10	....	2–1	832.5	0.38
	1021.32	....	2–2	832.5	1.12
C II	1036.330	$2p\ ^2P^0$–$2p^2\ ^2S$	$\frac{1}{2}$–$\frac{1}{2}$	0	0.24
	1037.017	....	$\frac{3}{2}$–$\frac{1}{2}$	64.0	0.48
S IV	1062.672	$3p\ ^2P^0$–$3p^2\ ^2D$	$\frac{1}{2}$–$\frac{3}{2}$	0	0.94
	1072.992	....	$\frac{3}{2}$–$\frac{5}{2}$	950.2	0.19
	1073.522	....	$\frac{3}{2}$–$\frac{3}{2}$	950.2	1.69
Cl II	1063.83	$3p^4\ ^3P$–$3p^5\ ^3P^0$	2–1	0	0.94
	1067.94	....	1–0	697	0.75
	1071.05	....	2–2	0	2.82
	1071.76	....	1–1	697	0.56
	1075.24	....	0–1	996	0.75
	1079.08	....	1–2	697	0.94
C II	1065.883	$2p^2\ ^2D$–$2p^3\ ^2P^0$	$\frac{5}{2}$, $\frac{3}{2}$–$\frac{3}{2}$	74932	0.87
	1066.121	....	$\frac{3}{2}$–$\frac{1}{2}$	74933.2	0.43
S III	1077.835	$3p^2\ ^1D$–$3p^3\ ^1D^0$	2–2	11320	3.50
N II	1083.977	$2p^2\ ^3P$–$2p^3\ ^3D^0$	0–1	0	0.67
	1084.568	....	1–1, 2	49.1	2.01
	1085.536	....	2–1, 2	131.3	0.54
	1085.699	....	2–3	131.3	2.82

TABLE 3—*Continued*

Ion	Wavelength (Å)	Multiplet	J_l–J_u	Excitation (cm^{-1})	gf
Si III	1108.35	$3p\ ^3P^0$–$3d\ ^3D$	0–1	52630	0.40
	1109.95	....	1–1, 2	52758	1.20
	1113.20	....	2–1, 2, 3	53019	2.00
Fe III	1122.526	$a\ ^5D$–$z\ ^5P^0$	4–3	0	4.50
	1124.883	....	3–2	436.2	2.32
	1126.72	....	2–1	738.9	0.88
	1128.02	....	3–3	436.2	1.17
	1128.72	....	2–2	738.9	1.46
	1129.19	....	1–1	932.4	1.12
	1130.404	....	0–1	1027.3	0.50
	1131.194	....	1–2	932.4	0.38
	1131.914	....	2–3	738.9	0.17
C III	1175.7	$2p\ ^3P^0$–$2p^2\ ^3P$	6 lines	52350	2.97
S III	1190.17	$3p^2\ ^3P$–$3p^3\ ^3D^0$	0–1	0	0.42
	1194.02	....	1–2	297.2	0.94
	1194.40	....	1–1	297.2	0.32
	1200.97	....	2–3	832.5	1.77
	1201.71	....	2–2	832.5	0.32
	1202.10	....	2–1	832.5	0.021
Si III	1206.52	$3s^2\ ^1S$–$3p\ ^1P^0$	0–1	0	1.68
C III	1247.368	$2p\ ^1P^0$–$2p^2\ ^1S$	1–0	102351.4	0.27
S II	1250.50	$3p^3\ ^4S^0$–$3p^4\ ^4P$	$\frac{3}{2}$–$\frac{1}{2}$	0	0.45
	1253.79	....	$\frac{3}{2}$–$\frac{3}{2}$	0	0.89
	1254.53	....	$\frac{3}{2}$–$\frac{5}{2}$	0	1.34
Si III	1294.55	$3p\ ^3P^0$–$3p^2\ ^3P$	1–2	52758	0.50
	1296.72	....	0–1	52630	0.40
	1298.90	....	{2–2, 1–1}	52900	1.80
	1301.12	....	1–0	52758	0.40
	1303.30	....	2–1	53019	0.50
C II	1323.916	$2p^2\ ^2D$–$2p^3\ ^2D^0$	4 lines	74932	1.7
C II	1334.515	$2p\ ^2P^0$–$2p^2\ ^2D$	$\frac{1}{2}$–$\frac{3}{2}$	0	0.92
	1335.648	....	$\frac{3}{2}$–$\frac{3}{2}$, $\frac{5}{2}$	64.0	1.84
Si IV	1393.73	$3s\ ^2S$–$3p\ ^2P^0$	$\frac{1}{2}$–$\frac{3}{2}$	0	1.31
	1402.73	....	$\frac{1}{2}$–$\frac{1}{2}$	0	0.65
C IV	1548.195	$2s\ ^2S$–$2p\ ^2P^0$	$\frac{1}{2}$–$\frac{3}{2}$	0	0.37
	1550.768	....	$\frac{1}{2}$–$\frac{1}{2}$	0	0.19
C II	1760.40	$2p^2\ ^2D$–$3p\ ^2P^0$	$\frac{5}{2}$, $\frac{3}{2}$–$\frac{3}{2}$	74932	3.9
	1760.81	....	$\frac{3}{2}$–$\frac{1}{2}$	74933.2	2.0
Al III	1854.722	$3s\ ^2S$–$3p\ ^2P^0$	$\frac{1}{2}$–$\frac{3}{2}$	0	1.31
	1862.782	....	$\frac{1}{2}$–$\frac{1}{2}$	0	0.65
Fe III	1895.456	$a\ ^7S$–$z\ ^7P^0$	3–4	30088.84	10.8
	1914.056	....	3–3	30088.84	8.4
	1926.304	....	3–2	30088.84	6.0
C III	2297.60	$2p\ ^1P^0$–$2p^2\ ^1D$	1–2	102351.4	0.72

Atomic Energy Levels (Moore, 1949). The gf-values for hydrogen were taken from the calculations of Underhill and Waddell (1959). All the rest were obtained from Gaustad and Spitzer (1961), Varsavsky (1961), or Allen (1963). Only a few of these were based on reliable measurements or calculations; most were estimated using sum rules. The partition functions of the ions were represented by the ground-state statistical weights.

III. COMPUTATION PROCEDURE

The effect of the blended line profiles on the emergent flux was computed using a program kindly provided by Miss Underhill (1962*b*). A set of 660 wavelengths was chosen from 911.6 to 3000 Å, spaced according to the irregularities of the spectrum. At each wavelength and at each of forty-four levels specified by a particular temperature and electron pressure, the absorption of each line was calculated from the abundance of the element, the populations of the ion state and the lower level, the *gf*-value, the broadening mechanism, and the distance from the line center. These opacities were summed over all contributing lines and added to the bound-free absorption of H I, H^-, He I, and He II and the scattering of the free electrons. An iterative procedure, described by Underhill (1960), was used to solve the transfer equation for the source function $S_\nu(t_\nu)$ at each wavelength and each depth t_ν. The monochromatic flux at the surface $F_\nu(0)$ was then found from the integral

$$F_\nu(0) = 2\int_0^\infty S_\nu(x)E_2(x)\,dx\,,$$

where $E_2(x)$ is the second exponential integral. No attempt was made to iterate the model to obtain flux constancy throughout the atmosphere.

The broadening in the Lyman lines of hydrogen was represented by a simplified expression derived by Griem (1960), which included the effects of ions by the quasi-static approximation and the effects of electrons by the impact approximation. Griem (1962) has since added two corrections at large separations from the line center—one a reduction because only the electron interaction times of shortest duration are significant, and the other an increase, even further in the wings, for the quasi-static effects of electrons. For all other lines a Doppler-damping profile was assumed with a damping constant ten times the classical value, so that

$$\delta = \frac{10}{4\pi}\frac{8\pi^2 e^2}{3m_e c}\frac{1}{\lambda_0^2} = \frac{0.177}{\lambda_0^2}\ \mathrm{sec}^{-1}\,.$$

Rudkjobing (1949) has noted that in B stars, where most of the hydrogen is ionized, electrons make the principal contribution to the damping, but no numbers are available for ultraviolet lines. The value adopted here is based on solar observations (Minnaert and Mulders 1931), although in the Sun the broadening is probably by the van der Waals forces of hydrogen atoms.

IV. DISCUSSION OF THE RESULTS

For the B2 V model, Figures 1–5 show the blended profiles of all the lines in Table 2 except Fe III at 1895 Å, which was left out to save space. It is only slightly wider than the other two in the multiplet in Figure 5. The continuum flux, without any lines, is represented by the nearly horizontal lines across the upper parts of the figures. The ordinates give fluxes F_ν in units of erg sec^{-1} cm^{-2} cps^{-1}, but a factor π is necessary to obtain the power radiated by the surface, so that

$$\sigma T_e^4 = \pi\int_0^\infty F_{\nu c}d\nu\,.$$

Several important features of the predicted spectrum are immediately apparent. The overlapping lines absorb one-half or more of the continuum flux over wide bands, particularly between 911.6 and 1040 Å, between 1060 and 1090 Å, and between 1165 and 1220 Å. The cores of all lines are very deep, a consequence of the low continuous opacity at these wavelengths and the low temperatures in the outermost layers. This characteristic is especially noticeable shortward of 1500 Å where the lines are on the steep side of the Planck function so that there is relatively little emission in the cores. The central depths are independent of abundance and *gf*-values but do decrease toward shorter wavelengths. This is consistent with the expression for the residual intensity at the core of a line given by Aller (1953) for an approximate solution of the transfer equation when the line is formed by pure absorption. Any scattering in the formation process would only

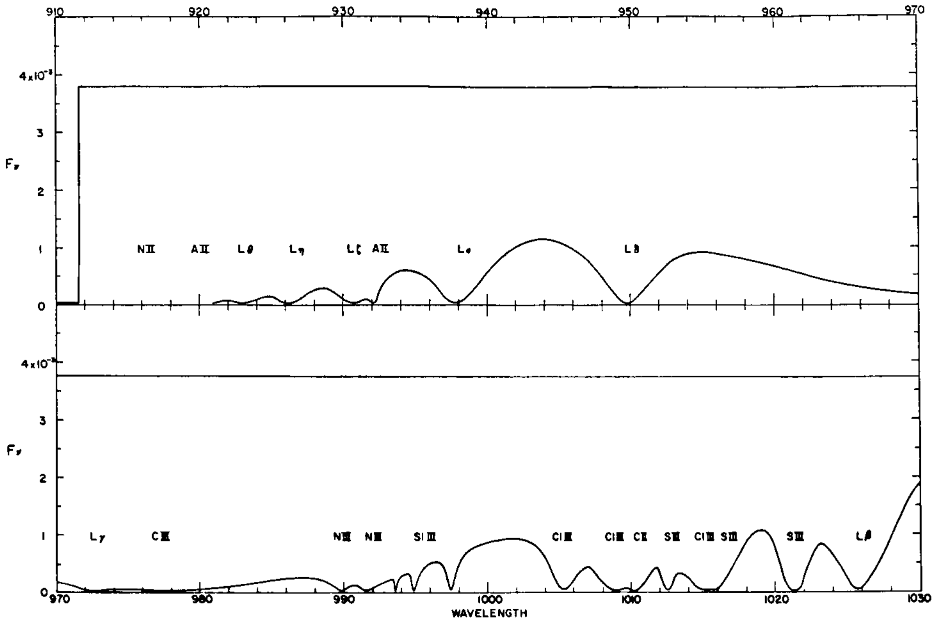

FIG. 1.—The spectrum of a B2 V star from 910 to 1030 Å

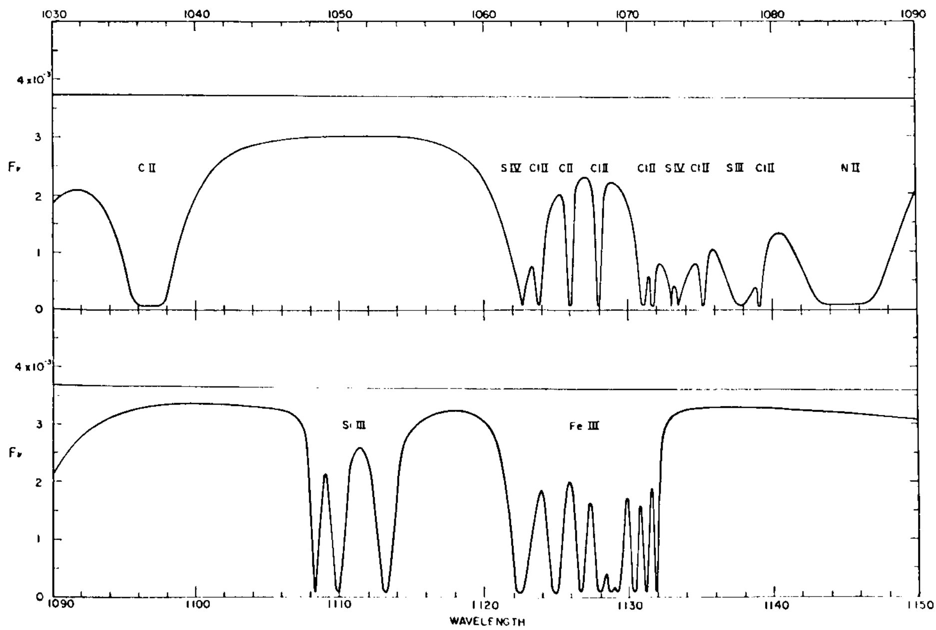

FIG. 2.—The spectrum of a B2 V star from 1030 to 1150 Å

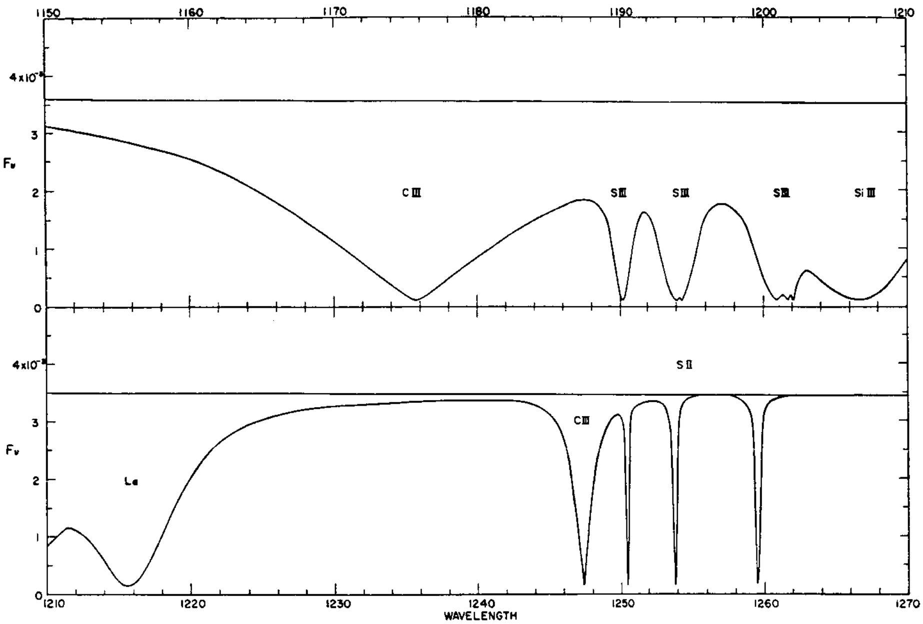

FIG. 3.—The spectrum of a B2 V star from 1150 to 1270 Å

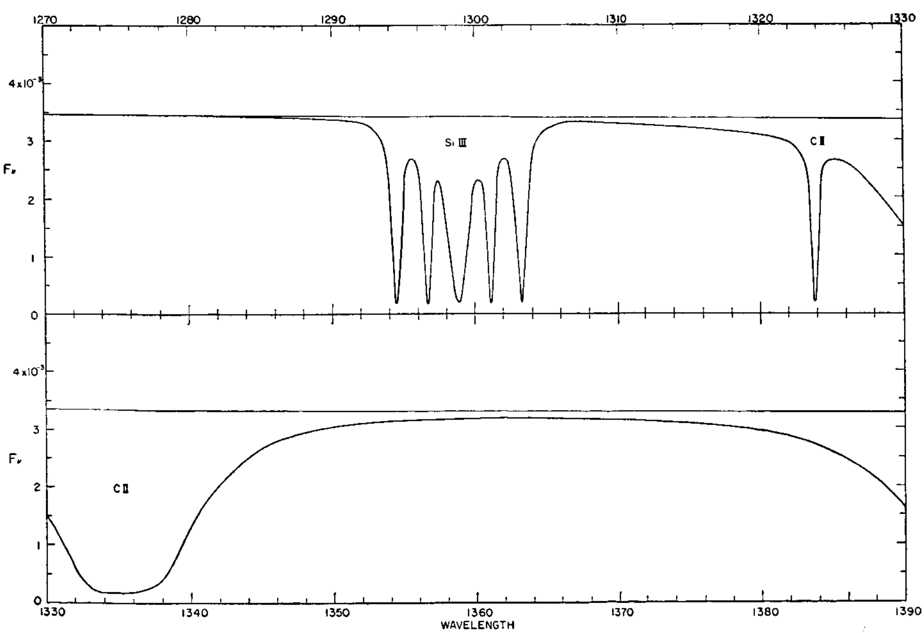

FIG. 4.—The spectrum of a B2 V star from 1270 to 1390 Å

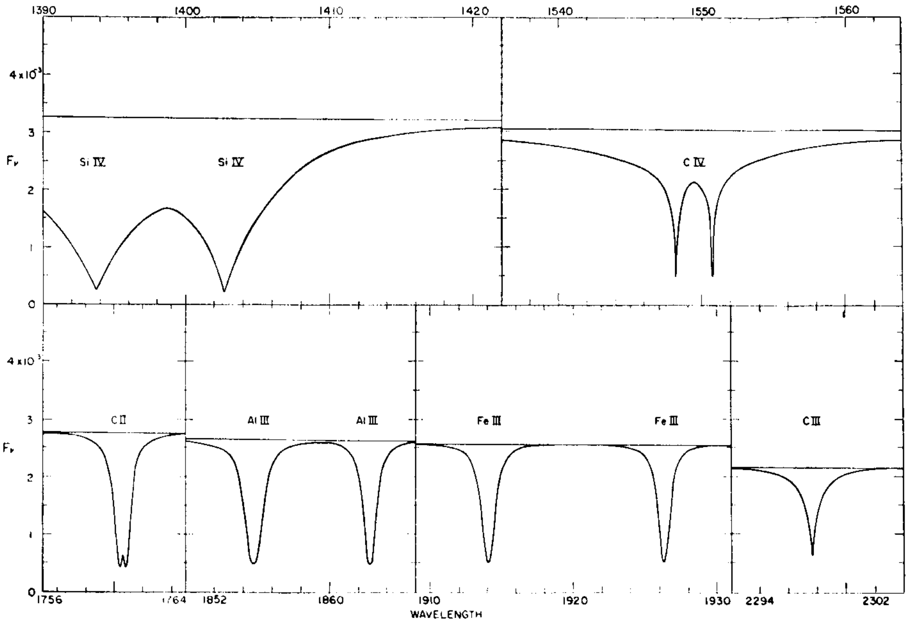

Fig. 5.—The spectrum of a B2 V star from 1390 to 3000 Å

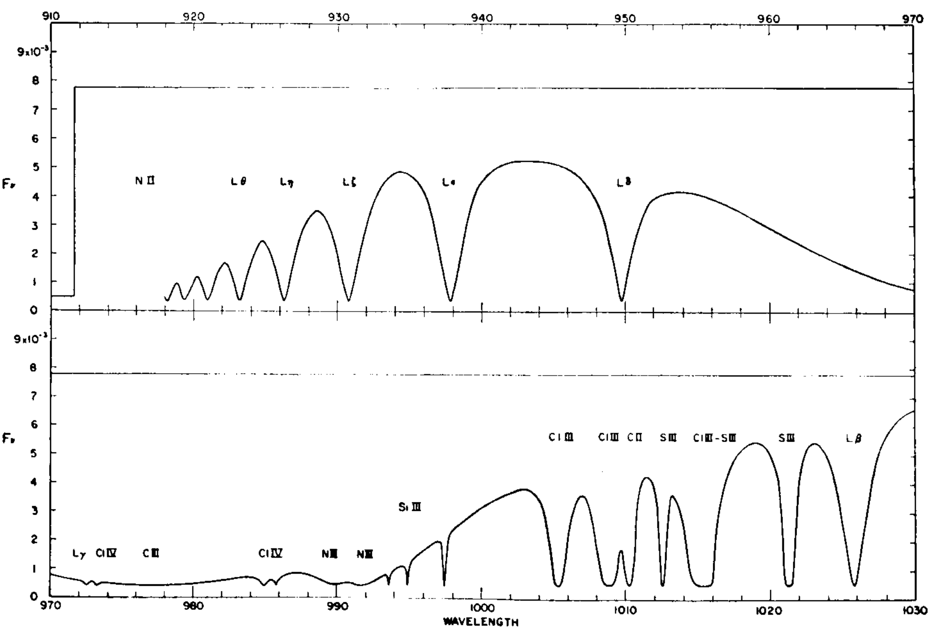

Fig. 6.—The spectrum of a B0 V star from 910 to 1030 Å

deepen a line further. Calculation of a few lines on the assumption that they were formed by pure scattering confirmed this and showed 10 per cent or less decrease in width. The precise values of the central intensities are not reliable, because at these points the opacity is so high that optical depth unity is reached in one or two integration steps. However for all fluxes greater than 0.3×10^{-3} there are sufficient steps.

Most of the lines shown are broader than any in the visual spectrum of a B star, where even the Balmer lines are only a few angstroms wide. These ultraviolet lines resemble more the H- and K-calcium lines in the Sun. The converging Lyman series dominates the spectrum shortward of 955 Å, and the equivalent widths increase a little with increasing quantum number. Longward of this wavelength the outstanding features are C III at 977.0 Å, N II at 1085 Å, C II at 1175.7 Å, Si III at 1206.5 Å, C II at 1335 Å, and Si IV at 1400 Å. All these lines are wider than 10 Å and the wings of the 977.0 line can be detected beyond 1050 Å. Below 1225 Å the continuum is seen in just a few narrow bands around 1050, 1100, 1118, and 1140 Å.

In Figure 6 only the first 120 Å of the B0 V spectrum longward of the Lyman limit is reproduced, but this is sufficient to show the general characteristics. Here the Lyman series and most of the other lines are somewhat weaker than in the B0 V model so that more of the continuum is visible. However, a few lines in the higher ionization states are stronger, such as N III at 990, S IV at 1062.7 and 1073, Si IV at 1400, and C IV at 1550. The central intensities are a little higher. The He II line at 1640 Å, corresponding to Hα, is only 1 Å wide.

TABLE 4

LINE BLANKETING

Spectrum	T_e (°K)	F_c (erg cm^{-2} sec^{-1})	$\Delta F/F_c$
B2 V......	25673	7.84×10^{12}	0.31
B0 V......	31023	1.67×10^{13}	0.20

The area between the two curves in the figures corresponds to the total flux absorbed by the lines. This correction of the calculated flux already has been applied in regions around 1314 and 1427 Å (Morton 1964) to reduce the discrepancies between the models and the ultraviolet measures of B stars with broad-band detectors in rockets. For the two models, Table 4 lists the effective temperature T_e, the integrated continuum flux F_c, and the fraction $\Delta F/F_c$ redistributed to other wavelengths by the lines. There is no major contribution to the line absorption longward of 3000 Å; the Balmer blanketing amounts to only 0.003 at B2 V and to one-half that at B0 V. In these calculations some strong lines may have been omitted; furthermore, Miss Underhill (1963) has suggested that many weak lines could contribute significantly in some parts of the spectrum. Therefore the values for $\Delta F/F_c$ quoted here must be considered as lower limits. Strömgren (1964) studied the effects of the ultraviolet lines in an approximate way by constructing models with the Lyman limit shifted 200 Å to longer wavelengths. For the same visual spectrum the effective temperature was lowered by about 2 per cent. Thus the blanketing should have little effect on the bolometric corrections.

V. SUMMARY

The ultraviolet spectra of two early B stars have been calculated from 911.6 to 3000 Å including the blended profiles of all absorption lines expected to be wider than 2 Å. The overlapping wings absorb a large fraction of the continuum flux shortward of 1410 Å producing a significant blanketing effect. At $T_e = 25673$, this amounts to 0.31 of the flux, while at $T_e = 31023$ it is 0.20. The former value is comparable with the fraction

0.29 found by Gaustad and Spitzer for an isothermal atmosphere at 20000° K. These results demonstrate the importance of including the opacities of the strong ultraviolet metal lines as well as the Lyman series when constructing a B-star atmosphere.

The details of the spectra shown here are necessarily questionable because the effects of the lines on the flux constancy have not been determined. Furthermore, since the temperatures in the outermost layers probably are too high and since these layers are not covered by enough integration steps, the cores of the lines are the last reliable. Additional uncertainties arise from the f-values and the damping constants used.

More accurate theoretical treatments can lead to improvements in all these areas, but other problems will be resolved only with measurements from above the atmosphere. They are necessary for information about abundances, chromospheric effects, the ratio of scattering to absorption in the line, and the contribution of weaker lines to the blanketing. Of course these observations may add many other features which have not been anticipated. However, this preliminary survey of the ultraviolet spectrum should be useful in planning instrumentation and observing programs for astronomical rockets and satellites. Gaustad and Spitzer already have noted that the possible interstellar line of molecular hydrogen at 1049 Å coincides with a continuum window in the B2 V star. The results for the hotter models suggest that more spectral regions clear of stellar lines will be found in earlier types.

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